arXiv:astro-ph/9911130v3 15 Sep 2000

Evolution of Neutron-Star, Carbon-Oxygen White-Dwarf Binaries

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ABSTRACT

At least one, but more likely two or more, eccentric neutron-star, carbon-oxygen white-dwarf binaries with an unrecycled pulsar have been observed. According to the standard scenario for evolving neutron stars which are recycled in common envelope evolution we expect to observe $\gtrsim 50$ such circular neutron star-carbon oxygen white dwarf binaries, since their formation rate is roughly equal to that of the eccentric binaries and the time over which they can be observed is two orders of magnitude longer, as we shall outline. We observe at most one or two such circular binaries and from that we conclude that the standard scenario must be revised.

Introducing hypercritical accretion into common envelope evolution removes the discrepancy by converting the neutron star into a black hole which does not emit radio waves, and therefore would not be observed.

Subject headings: binaries: close – stars: neutron – white dwarfs – stars: evolution – stars: statistics

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1. Introduction

We consider the evolution of neutron-star, carbon-oxygen white-dwarf binaries using both the Bethe & Brown (1998) schematic analytic evolutions and the Portegies Zwart & Yungelson (1998) numerical population syntheses.

The scenario in which the circular neutron-star, carbon-oxygen white-dwarf binaries (which we denote as $(ns, co)_{\mathcal{C}}$ hereafter) have gone through common envelope evolution is considered. In conventional common envelope evolution for the circular binaries it is easy to see that the observed ratio of these to eccentric binaries (hereafter $(ns, co)_{\mathcal{E}}$) should be ~ 50 because: (i) The formation rate of the two types of binaries is, within a factor 2, the same. (ii) The magnetic fields in the circular binaries will be brought down by a factor of ~ 100 by He accretion in the neutron-star, He-star phase following common envelope evolution just as the inferred pulsar magnetic field strengths in the double neutron star binaries are brought down (Brown 1995). In the eccentric binaries the neutron star is formed last, after the white dwarf, so there is nothing to circularize its orbit. More important, its magnetic field will behave like that of a single star and will not be brought down from the $B \sim 10^{12}$ G with which it is born. (At least empirically, neutron star magnetic fields are brought down only in binaries, by accreting matter from the companion star, Taam & Van den Heuvel 1986, although Wijers 1997 shows the situation to be more complex.) Neutron stars with higher magnetic fields can be observed only for shorter times, because of more rapid spin down from magnetic dipole radiation. The time of possible observation goes inversely with the magnetic field B. We use the observability premium

$$\Pi = 10^{12}\mathrm{G}/B \tag{1}$$

(Wettig & Brown 1996) which gives the relative time a neutron star can be observed. Given our above point (ii), the circular binaries should have an observability premium $\Pi \sim 100$ as compared with $\Pi \sim 1$ for the higher magnetic field neutron star in an eccentric orbit. Correcting for the factor 2 higher formation rate of the eccentric binaries (point (i) above) this predicts the factor ~ 50 ratio of circular to eccentric binaries.

In our paper we cite one firm eccentric neutron-star, carbon-oxygen white-dwarf binary $(ns, co)_{\mathcal{E}}$ B2303+46 and argue for a recently observed second one, J1141−65. Portegies Zwart & Yungelson (1999) suggest PSR 1820−11 may also be in this class, but cannot exclude the possibility that the neutron star companion is a main sequence star (Phinney & Verbunt 1991). This would imply that $\gtrsim 100$ such binaries with circular orbits should be observed. But, in fact, only one [2] B0655+64 is observed if we accept the developing

[2] We have a special scenario for evolving it; see section 3.2.

concensus (Section 2.3) that those observed $(ns, co)_{\mathcal{C}}$ are evolved with avoidence of common envelope evolution. We are thus confronted by a big discrepancy, for which we suggest a solution.

In order to understand our solution, we need to review three past works. In the earlier literature the observed circular $(ns, co)_{\mathcal{C}}$ were evolved through common envelope, e.g., see Van den Heuvel (1994) and Phinney & Kulkarni (1994). Accretion from the evolving giant progenitor of the white dwarf was neglected, since it was thought that the accretion would be held to the Eddington rate of $\dot{M}_{\rm Edd} \sim 1.5 \times 10^{-8}\ M_\odot\ {\rm yr}^{-1}$, and in the ~ 1 year long common envelope evolution a negligible amount of matter would be accreted. We term this the standard scenario. However, Chevalier (1993) showed that once $\dot{M}$ exceeded $\sim 10^4 \dot{M}_{\rm Edd}$, it was no longer held up by the radiative pressure due to the X rays from the neutron star, but that it swept them inwards in an adiabatic inflow. Bethe & Brown (1998) employed this hypercritical accretion in their evolution of double neutron star (ns, ns) and neutron-star, low-mass black-hole binaries $(ns, lmbh)$ and we shall use the same techniques in binary evolution here. In particular, these authors found that including hypercritical accretion in the standard scenario for double-neutron star binary evolution, the first born neutron star went into a low-mass black hole. To avoid the neutron star going through the companion's envelope, a new scenario was introduced beginning with a double He star binary. It gives about the right number of double neutron star binaries.

A new development has been that most of the circular $(ns, co)_{\mathcal{C}}$ are currently evolved with avoidance of common envelope evolution. In Section 2.3, we shall summarize this work, carried out independently by King & Ritter (1999) and Tauris, Van den Heuvel & Savonije (2000). If we accept the new scenario, at most one or two circular $(ns, co)_{\mathcal{C}}$ that went through common envelope evolution have been observed. Yet, in the standard scenario at least ~ 50 of them should be seen.

In this paper we find that in (ns, co) which do go through common envelope evolution, the neutron star goes into a black hole. The (ns, co) binaries observed to date have been identified through radio emission from the neutron star. Thus, binaries containing a low-mass black hole would not have been seen. We discuss masses for which neutron stars go into black holes.

Although the main point of our paper relies only on relative formation rates, we shall show in Appendix A that the Bethe & Brown (1998) schematic, analytic analysis agrees well with the detailed numerical population synthesis of Portegies Zwart, once both are normalized to the same supernova rate.

2. The Problem

2.1. Standard Scenario vs Observations

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Portegies Zwart & Yungelson (1998), in a very careful population synthesis, have calculated the expected number of newly created binaries of compact stars (neutron stars or black holes) and white dwarfs. Among the latter, they distinguish between those consisting of helium and those consisting of carbon-oxygen (denoted *co*). To make an eccentric binary containing a neutron star, the supernova must occur after the carbon-oxygen star has formed (Portegies Zwart & Yungelson 1999). To make a circular binary containing a neutron star it is necessary that its companion be close so that at some earlier stage in evolution (but after formation of the neutron star) there was mass transfer or strong tidal interaction, which requires the companion to (nearly) fill its Roche Lobe.

Since the Bethe & Brown (1998) schematic calculations did not include mass exchange, which is very important in evolving $(ns, co)_{\mathcal{E}}$ binaries, we need the more complete calculations of Portegies Zwart & Yungelson, which are listed in Table 1. We discuss this in more detail later. These do not include hypercritical accretion; i.e., they follow the standard scenario. In this case the formation ratio of $(ns, co)_{\mathcal{C}}$ to $(ns, co)_{\mathcal{E}}$ is $17.7/32.1 = 0.55$. We now make the case that if the $(ns, co)_{\mathcal{C}}$ were to be formed through common envelope evolution (Phinney & Kulkarni 1994; Van den Heuvel 1994) in the standard scenario, their pulsar magnetic fields would be brought down to $B \sim 10^{10}$ G because of the similarity to binary neutron star systems in which this occurs. In detail, this results from helium accretion during the neutron-star, He-star stage which precedes the final binary (Brown 1995, Wettig & Brown 1996).

Detailed calculation of Iben & Tutukov (1993) for original donor masses $4-6\ M_\odot$ of the white dwarf progenitor show that following common envelope evolution the remnant stars fill their Roche lobes and continue to transfer mass to their companion neutron star. These remnants consist of a degenerate carbon-oxygen core and an evolving envelope undergoing helium shell burning. The mass transfer to the neutron star is at a rate $\dot{M} < 10^4 \dot{M}_{\rm Edd}$, the lower limit for hypercritical accretion, so it is limited by Eddington. Van den Heuvel (1994) estimates that the neutron star accretes about $0.045\ M_\odot$ and $0.024\ M_\odot$ in the case of the ZAMS $5\ M_\odot$ and $6\ M_\odot$ stars, and $0.014\ M_\odot$ for a $4\ M_\odot$ star, where these ZAMS masses refer to the progenitors of the white dwarfs. The accretion here is of the same order, roughly double,[3] the wind accretion used by Wettig & Brown (1996) in the evolution of

[3] The He burning time to be used for the progenitor of the white dwarf is $\sim 10^6$ years, whereas for the

the relativistic binary pulsars B1534+12 and B1913+16. There the magnetic fields were brought down by a factor ~ 100 from $B \sim 10^{12}$ G to $\sim 10^{10}$ G, increasing the observability premium Π by a factor of ~ 100. Thus, the scenario in which the $(ns, co)_{\mathcal{C}}$ are produced through common envelope evolution without hypercritical accretion should furnish them with $\Pi \sim 100$, by helium accretion following the common envelope. Although the detailed description may not be correct, the similarity of evolution of $(ns, co)_{\mathcal{C}}$ with that of binary neutron stars in the older works (Phinney & Kulkarni 1994; Van den Heuvel 1994) should furnish these with about the same Π.

There is one confirmed $(ns, co)_{\mathcal{E}}$, namely B2303+46, see Table 2, so there should be about 50 circular ones which went through common envelope evolution. Indeed, several circular ones have been observed (see Table 2), and one or two of these may have gone through common envelope evolution. Thus we have a big discrepancy between the standard scenario and the observations. In the next section, we discuss the possibility of PSR J1141−6545 being $(ns, co)_{\mathcal{E}}$, which enhances the discrepancy.

2.2. Is PSR J1141−6545 $(ns, co)_{\mathcal{E}}$?

Not only is the eccentric B2303+46 quite certain, but a relativistic counterpart, PSR J1141−6545 has recently been observed (Kaspi et al. 2000), in an eccentric orbit. The inferred magnetic dipole strength is 1.3×10^{12} G, and the total mass is $2.300 \pm 0.012\ M_{\odot}$. Kaspi et al. argue that the companion of the neutron star can only be a white dwarf, or neutron star. With a total mass of 2.3 $M_{\odot}$, if J1141−65 were to contain two neutron stars, each would have to have a mass of $\sim 1.15\ M_{\odot}$, well below the 19 accurately measured neutron star masses, see Fig. 1 (Thorsett & Chakrabarty 1999).

We can understand the absence of binary neutron stars with masses below $\sim 1.3\ M_{\odot}$, although neutron stars of this mass are expected to result from the relatively copious main sequence stars of ZAMS mass $\sim 10 - 13\ M_{\odot}$ from the argument of Brown (1997). The He stars in the progenitor He-star, pulsar binary of mass $\lesssim 4\ M_{\odot}$ (Habets 1986) expand substantially during He shell burning. Accretion onto the nearby pulsar sends it into a black hole. Indeed, with inclusion of mass loss by helium wind, He stars of masses up to 6 or 7 $M_{\odot}$ expand in this stage (Woosley, Langer & Weaver 1995). Fryer & Kalogera (1997) find that special kick velocities need to be selected in order to avoid the evolution of PSR 1913+16 and PSR 1534+12 from going into a black hole by reverse Case C mass

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relativistic binary pulsars the average time of 5×10^5 years is more appropriate, so one would expect a factor ~ 2 greater accretion.

transfer (mass transfer from the evolving He star companion onto the pulsar in the He-star, neutron-star stage which precedes that of the binary of compact objects).

Our above argument says that the first neutron star formed in these would be sent into a black hole when its companion He star evolved and poured mass on it. Therefore, we believe the companion in J1141−65 must be a white dwarf. Earlier Tauris & Sennels (2000) developed the case that J1141−65 was an eccentric neutron-star, white-dwarf binary. Given the high magnetic field of J1141−65 (1.3×10^{12} G) with low observability premium of 0.77, this would increase the predicted observed number of circular $(ns, co)_{\mathcal{C}}$ which had gone through common envelope evolution to ~ 130 in the standard scenario.

2.3. Evolution of Neutron-Star, Carbon-Oxygen White-Dwarf Binaries with Avoidance of Common Envelope Evolution

Our discussion of the common envelope evolution in the last section applied to convective donors. In case the donor is radiative or semiconvective, common envelope evolution can be avoided. Starting from the work of Savonije (1983), Van den Heuvel (1995) proposed that most low mass X-ray binaries would evolve through a Her X-1 type scenario, where the radiative donor, more massive than the neutron star, poured matter onto its accretion disk at a super Eddington rate, during which time almost all of the matter was flung off. This involved Roche Lobe overflow. Although Van den Heuvel limited the ZAMS mass of the radiative donor to 2.25 $M_\odot$ in order to evolve helium white-dwarf, neutron star binaries, his scenario has been extended to higher ZAMS mass donors in order to evolve the carbon-oxygen white-dwarf, neutron star binaries. The advection dominated inflow-outflow solutions (ADIOS) of Blandford & Begelman (1999) suggest that the binding energy released at the neutron star can carry away mass, angular momentum and energy from the gas accreting onto the accretion disk provided the latter does not cool too much. In this way the binding energy of gas at the neutron star can carry off $\sim 10^3$ grams of gas at the accretion disk for each gram accreting onto the neutron star. King & Begelman (1999) suggest that such radiatively-driven outflows allow the binary to avoid common envelope evolution.

As noted above, for helium white dwarf companions, Van den Heuvel (1995) had suggested Cyg X-2 as an example following the Her X-1 scenario. King & Ritter (1999) calculated the evolution of Cyg X-2 in the ADIOS scenario in detail. These authors also evolved the $(ns, co)_{\mathcal{C}}$ binaries in this way, using donor stars of ZAMS masses $4 - 7$ $M_\odot$. Tauris, Van den Heuvel, & Savonije (1999) have carried out similar calculations, with stable mass transfer. These authors find that even for extremely high mass-transfer rates, up to

$\dot{M} \sim 10^4 \dot{M}_{\rm Edd}$, the system will be able to avoid a common envelope and spiral-in evolution.

Tauris, van den Heuvel & Savonije 2000 evolve J1453−58, J1435−60 and J1756−5322, the three lowest entries in our Table 2, through common envelope. We obtained the eccentricities and $\dot{P}$'s for the first two of these (Fernando Camilo, private communication). The binary J1453−58, quite similar to J0621+1002 has a substantial eccentricity and clearly should be evolved with a convective donor as Tauris et al did for J0621+1002. The spin periods of J1435−60 and J1756−5322 are short, indicating greater recycling than the other listed pulsars. It would seem difficult to get the inferred magnetic field down to the 4.7×10^8 G of J1435−60 by the Iben & Tutukov or Wettig & Brown accretion scenarios following common envelope evolution as discussed in Section 2.1. If, however, one does believe that J1435−60 and J1756−5322 have gone through common envelope, the discrepancy between predicted and observed circular binaries in the standard scenario is only slightly relieved.

2.4. Are There Observational Selection Effects ?

In Table 3 we have tabulated $S_{400} \times d^2$ in order to see whether the normalized intensity gives strong selection effects. Note that the 35.95 for B2303+46 is not so different from the 43.56 and 203.35 for B1534+12 and B1913+16, respectively. For the circular binaries $(ns, co)_{\mathcal{C}}$ the intensities are less, but their empirical Observability Premium Π is much larger. There may be other observational selection effects, but, we believe that there are no observational selection effects strong enough to compensate for the factor 100 discrepancy between the observed population and the one expected from the standard model. So the problem remains the same.

3. The Answer

3.1. Black Hole Formation in Common Envelope Evolution

We believe the answer to the missing binaries is that the neutron star goes into a black hole in common envelope evolution, as we now describe. We label the mass of the neutron star as M_A and that of the giant progenitor of the white dwarf as M_B. Following Bethe & Brown (1998) we choose as variables the neutron star mass M_A and $Y \equiv M_B/a$, where a is the orbital radius. From their eq. (5.12) we find

$$\frac{M_{A,f}}{M_{A,i}} = \left(\frac{Y_f}{Y_i}\right)^{c_d - 1} \qquad (2)$$

where c_d is the drag coefficient. From Shima et al. (1985) we take

$$c_d = 6. \tag{3}$$

We furnish the energy to remove the hydrogen envelope of the giant B (multiplied by α_{ce}^{-1}, where α_{ce} is the efficiency of coupling of the orbital motion of the neutron star to the envelope of B) by the drop in orbital energy of the neutron star; i.e.,

$$\frac{0.6\ GM_{B,i}Y_i}{\alpha_{ce}} = \frac{1}{2}GM_{A,i}Y_i\left(\frac{Y_f}{Y_i}\right)^{6/5}. \tag{4}$$

Here the $0.6GM_{B,i}Y_i$ is just the binding energy of the initial giant envelope, found by Applegate (1997) to be $0.6GM_{B,i}^2a_i^{-1}$, and the right hand side of the equation is the final gravitational binding energy $\frac{1}{2}GM_{A,f}M_{B,f}a_f^{-1}$ in our variables. Using eqs. (2) and (3) in eq. (4) one finds

$$\frac{M_{A,f}}{M_{A,i}} = \left(\frac{1.2M_{B,i}}{\alpha_{ce}M_{A,i}}\right)^{1/c_d}. \tag{5}$$

For the sake of argument, we take the possible range of initial neutron star mass to be $1.2 - 1.5\ M_\odot$ (the upper bound is the Brown & Bethe (1994) mass at which a neutron star goes into a low-mass black hole), and the main sequence progenitor masses of the carbon-oxygen white dwarf to be $M_{B,i} = 2.25 - 10\ M_\odot$. As we show in Appendix C, in the Bethe & Brown (1998) schematic model, mass transfer was assumed to take place when the evolving giant reached the neutron star, whereas more correctly it begins when the envelope of the giant comes to its Roche Lobe. For the masses we employ, main sequence progenitors of the carbon-oxygen white dwarf of $2.25 - 10\ M_\odot$, the fractional Roche Lobe radius is

$$r_L \sim 0.5. \tag{6}$$

The binding energy of the progenitor giant at its Roche Lobe is, thus, double what it would be at a_i, the separation of giant and neutron star. Therefore, a Bethe & Brown $\alpha_{ce} = 0.5$ corresponds to a true efficiency $\hat{\alpha}_{ce} \sim 1$, if the latter is defined as the value for which the envelope removal energy, at its Roche Lobe, is equal to the drop in neutron star orbital energy as it moves from a_i to a_f. If we take $\alpha_{ce} = 0.5$ in eq. (5) we find, given our assumed possible intervals

$$1.54\ M_\odot \lesssim M_{A,f} \lesssim 2.38\ M_\odot. \tag{7}$$

These are above the neutron star mass limit $1.5\ M_\odot$ (Brown & Bethe 1994) beyond which a neutron star goes into a low-mass black hole. Thus, all neutron stars with common envelope evolution in our scenario evolve into black holes. This solves the big discrepancy between the standard scenario and observation. The only remaining problem is the evolution of B0655+64, which survived the common envelope evolution, and we suggest a special scenario for it in the next section.

3.2. Is B0655+64 a problem?

Van den Heuvel & Taam (1984) were the first to notice that the $(ns, co)_{\mathcal{C}}$ system B0655+64 might have been formed in a similar way as the double neutron stars. The short period of 1.03 days, magnetic field $\sim 10^{10}$ G, and the high companion mass of $\sim 1\ M_\odot$ make this binary most similar to a binary neutron star, but with a carbon-oxygen white-dwarf companion, resulting from probable ZAMS masses $\sim 5-8\ M_\odot$. For a 1.4 $M_\odot$ neutron star with 1 $M_\odot$ white-dwarf companion $a_f = 5.7 R_\odot$.

The similarity of B0655+64 to the close neutron star binaries suggests the double helium star scenario (Brown 1995) to calculate the evolution. The ZAMS mass of the primary is chosen to be just above the limit for going into a neutron star, that of the secondary just below. For the double He star scenario the ZAMS masses of primary and secondary cannot be more than $\sim 5\%$ different. However, in this case the ratio q of masses is so close to unity that the secondary will not be rejuvenated (Braun & Langer 1995: If the core burning of hydrogen to helium in the companion star is nearly complete, the accreted matter would have to cross a molecular weight barrier in order to burn and if q is near unity there is not time enough to do so. Thus He cores of both stars will evolve as if the progenitors never had more than their initial ZAMS mass.)

What we have learned recently about effects of mass loss (Wellstein & Langer 1999) will change the Brown (1995) scenario in detail, but not in general concept. An $\sim 10\ M_\odot$ ZAMS star which loses mass in RLOF to a lower mass companion will burn helium as a lower-mass star due to subsequent mass loss by helium winds, roughly as an 8 $M_\odot$ star (Wellstein & Langer, in preparation). Thus, the primary must have ZAMS mass $\gtrsim 10\ M_\odot$ in this case in order to evolve into a neutron star following mass loss. Although the secondary will not be rejuvenated as mass is transferred to it, it will burn helium without helium wind loss because it is clothed with a hydrogen envelope. Thus, a secondary of ZAMS 8 $M_\odot$ will burn He roughly as the primary of 10 $M_\odot$ in the situation considered. Given these estimates, a primary of ZAMS mass $M \lesssim 10\ M_\odot$ will evolve into a white dwarf, whereas a secondary of mass $\gtrsim 8\ M_\odot$ will end up as a neutron star. Of course, the former must be more massive than the latter, but stars in this mass range are copious because this is the lowest mass range from which neutron stars can be evolved, so there will be many such cases.

This scenario might not be as special as outlined because the fate of stars of ZAMS mass $8-10\ M_\odot$, which do not form iron cores but do burn in quite different ways from more massive stars, is somewhat uncertain in the literature. Whereas it is generally thought that single stars in this range end up as neutron stars, it has also been suggested that some of them evolve as AGB stars ending in white dwarfs. In terms of these discussions it does not seem unlikely that with two stars in the binary of roughly the same mass, the first

to evolve will end up as a neutron star and the second as a white dwarf, especially if the matter transferred in RLOF cannot rejuvenate the companion.

Van den Heuvel & Taam (1984) evolved B0655+64 by common envelope evolution. In taking up the problem again, Tauris, Van den Heuvel & Savonije (2000) in agreement with King & Ritter (1999) find that B0655+64 cannot be satisfactorily evolved with their convective donor scenario. Tauris et al. suggest a spiral-in phase is the most plausible scenario for the formation of this system, but we find that the neutron star would go into a black hole in this scenario, unless the two progenitors burn He at the same time.

3.3. Neutron Star Masses

There is by no means agreement about maximum and minimum neutron star masses in the literature. The mass determination of Vela X-1 have been consistently higher than the Brown & Bethe 1.5 $M_\odot$ which is consistent with well measured neutron star masses in Fig. 1. In a recent careful study at ESO Barziv et al. (2000), as reported by Van Kerkwijk (2000), obtain

$$M_{NS} = 1.87^{+0.23}_{-0.17} \ \ M_\odot. \tag{8}$$

Even at 99% confidence level, $M_{NS} > 1.6\ M_\odot$. Taking the maximum mass to be 1.87 $M_\odot$ and $\alpha_{ce} = 0.5$, $(M_{NS})_{min} = 1.2\ M_\odot$ one finds from eq. (5) that the maximum carbon-oxygen white dwarf progenitor mass of $(ns, co)_{\mathcal{C}}$ is

$$(M_{B,i})_{max} = \alpha_{ce} \left(\frac{M_{A,f}}{M_{A,i}}\right)^6 \frac{M_{A,i}}{1.2} \approx 7.2\ M_\odot. \tag{9}$$

Although there is some uncertainty in the efficiency α_{ce}, the ratio $M_{B,i}/\ M_\odot$ is much more sensitive to $M_{A,f}$ because of the 6th power of the ratio in eq. (9).[4] But then one cannot explain why no $(ns, co)_{\mathcal{C}}$ (except B0655+64) which survived the common envelope evolution are seen, since this mass is high enough to give those of the white dwarf companions.

Distortion of the $\sim 20\ M_\odot$ B-star companion by the neutron star in Vela X-1 brings in large corrections (Zuiderwijk et al. 1977, van Paradijs et al. 1977a, van Paradijs et al. 1997b) making measurement of neutron star masses in high-mass X-ray binaries much more difficult than those with degenerate companions.

[4] With $M_{NS} = 1.5\ M_\odot$ we get $(M_{B,i})_{max} = 1.9\ M_\odot$ which is below the minimum M_B ($\sim 2.25\ M_\odot$) for forming a carbon-oxygen white dwarf, so no $(ns, co)_{\mathcal{C}}$ survive the common envelope evolution.

Given the $Y_e \simeq 0.43$ at the collapse of the core of a large star (Aufderheide et al. 1990) one finds the cold Chandrasekhar mass to be

$$M_{CS} = 5.76\ Y_e^2\ M_\odot \approx 1.06\ M_\odot \tag{10}$$

where Y_e is the ratio of the number of electrons to the number of nucleons. Thermal corrections increase this a bit, whereas lattice corrections on the electrons decrease it, so that when all is said and done, $M_{CS} \gtrsim 1.1\ M_\odot$ (Shapiro & Teukolsky 1983). The major dynamical correction to this is from fallback following the supernova explosion. We believe that fallback in supernova explosions will add at least $\sim 0.1\ M_\odot$ to the neutron star, since bifurcation of the matter going out and in happens at about 4000 km (Bethe & Brown 1995). Thus our lower limit of $\sim 1.2\ M_\odot$ is reasonable.

4. Discussion and Conclusions

At least one, but more likely two or more, $(ns, co)_\mathcal{E}$ binaries with an unrecycled pulsar have been observed. According to the standard scenario for evolving neutron stars which are recycled in a common envelope evolution we then expect to observe $\gtrsim 50$ $(ns, co)_\mathcal{C}$. We only observe B0655+64 (which we evolve in our double He-star way) and possibly one or two binaries that went through common envelope evolution and from that we conclude that the standard scenario must be revised. Introducing hypercritical accretion into common envelope evolution (Brown 1995; Bethe & Brown 1998) removes the discrepancy.

We believe that the evolution of the other $(ns, co)_\mathcal{C}$ binaries may originate from systems with a neutron star with a radiative or semi-convective companion. The accretion rate in these systems can be as high as $10^4 \dot{M}_{\rm Edd}$ but common envelope evolution is avoided. This possibility, however, does not affect our conclusion concerning hypercritical accretion.

It is difficult to see "fresh" (unrecycled) neutron stars in binaries because they don't shine for long. B2303+46 (Table 2) is the most firm example of a $(ns, co)_\mathcal{E}$ binary with a fresh neutron star. Although binaries where a "fresh" neutron star is accompanied by a black hole have similar birthrates ($\sim 10^{-4}$ yr^{-1} for both types; Bethe & Brown, 1999, and Portegies Zwart & Yungelson 1999) and lifetime, none are observed. In Appendix A we quote results of Ramachandran & Portegies Zwart (1998) that show there is an observational penalty which disfavors the observation of neutron stars with black holes as companions, because of the difficulty in identifying the pulsar due to the Doppler shift which smears out the signal in these short-period objects. Because of the longer orbital period and lower companion mass of $(ns, co)_\mathcal{E}$, such binaries are less severely plagued by this effect, although the recently discovered J1141−6545 is a relativistic binary with 5 hr

period. We therefore argue that it is not unreasonable that no $(lmbh, ns)$ binaries have yet been observed, but that they should be actively searched for since the probability of seeing them is not far down from that of seeing the $(ns, co)_{\mathcal{E}}$'s.

We are grateful to Fernando Camilo for information in Table 2. We would like to thank Brad Hansen, Marten van Kerkwijk, Ralph Wijers, Thomas Tauris and Lev Yungelson for useful discussions and advice, and thank Justin Holmer for providing us with the convective envelopes for stars in the giant phase. GEB would like to thank Brad Hansen also for correspondence which started him on this problem. GEB & CHL were supported by the U.S. Department of Energy under grant DE-FG02-88ER40388.

SPZ was supported by NASA through Hubble Fellowship grant HF-01112.01-98A by the Space Telescope Science Institute, which is operated by the Association of Universities for Research in Astronomy, Inc., for NASA under contract NAS 5-26555. SPZ is grateful for the hospitality of the State University of New York at Stony Brook.

A. Comparison of Population Syntheses

In this Appendix we first compare results that the Bethe & Brown (1998) schematic analytic evolution would have given without hypercritical accretion with the Portegies Zwart & Yungelson (1998, 1999) results of Table 1, which do not include hypercritical accretion. We can then illustrate how hypercritical accretion changes the results.

Without hypercritical accretion the $(lmbh, ns)$ binaries of Bethe & Brown would end up rather as neutron star binaries (ns, ns), giving a summed formation rate of 1.1×10^{-4} yr^{-1}, to compare with 1.1×10^{-4} yr^{-1} from the Portegies Zwart numerical driven population sythesis results presented in Table 1. This good agreement indicates that kicks that the neutron star receives in formation were implemented in the same way in the two syntheses. Introduction of hypercritical accretion leaves only those neutron stars which do not go through a common envelope; i.e., those in the double He star scenario of Brown (1995), with formation rate 10^{-5} yr^{-1}. This is much closer to the estimated empirical rate of 8×10^{-6} yr^{-1} of Van den Heuvel & Lorimer (1996) which equals the rate derived independently by Narayan et al. (1991) and Phinney (1991). Large poorly known factors are introduced in arriving at these "empirical" figures, so it is useful that our theoretical estimates end up close to them. In our theoretical estimates the possibility described earlier that the pulsar in the lower mass binary pulsars goes into a black hole in the He shell burning stage of the progenitor He-star, neutron-star binary (Brown 1997) was not taken into account and this process may change $\sim$ half of the remaining neutron star binaries in our evolution into

$(lmbh, ns)$ binaries.

Bethe & Brown (1998) had a numerical symmetry between high-mass binaries in which both massive stars go supernova and those in which the more massive one goes supernova and the other, below the assumed dividing mass of 10 $M_\odot$, did not; i.e., the number of binaries was equal in the two cases. Taking ZAMS mass progenitors of $2.3 - 10\ M_\odot$ for carbon-oxygen white dwarfs, we then find a rate of

$$R = 2 \times \frac{10 - 2.3}{10} \times 1.1 \times 10^{-4}\ \mathrm{yr}^{-1} = 16.9 \times 10^{-5}\ \mathrm{yr}^{-1} \tag{A1}$$

for the formation rate of $(ns, co)_{\mathcal{C}}$ binaries. The 1.1×10^{-4} yr^{-1} is taken from the last paragraph and applies here because of the numerical symmetry mentioned above. The factor 2 results because there is no final explosion of the white dwarf to disrupt the binary, as there was above in the formation of the neutron star. This rate R is to be compared with $(ns, co)_{\mathcal{C}} = 17.7 \times 10^{-5}$ yr^{-1} in Table 1. These $(ns, co)_{\mathcal{C}}$ binaries are just the ones in which the neutron star goes into a black hole in common envelope evolution, unless the masses of the two initial progenitors are so close that they burn He at the same time. Then a binary such as B0655+64 can result, since the two helium stars then go through a common envelope, rather than the neutron star and main sequence star.

It is of interest to compare the populations of (ns, ns) binaries with the $(ns, co)_{\mathcal{E}}$ binaries. We must rely on the Portegies Zwart result for the latter, which cannot be evolved without mass transfer, which is not included in the Bethe & Brown evolution. The (ns, ns) binaries involve common envelope evolution, whereas the $(ns, co)_{\mathcal{E}}$ do not. Thus, results for the rates should differ substantially in the standard scenario, which does not include hypercritical accretion, and our scenario which does. The ratio for (ns, ns) and $(ns, co)_{\mathcal{E}}$ from Table 1 are 10.6×10^{-5} yr^{-1} and 32.1×10^{-5} yr^{-1}. The (ns, ns) are recycled in the He-star, pulsar stage by the He wind, giving an observability premium of $\Pi \sim 100$ (Brown 1995). The pulsar in the $(ns, co)_{\mathcal{E}}$ is not recycled. Thus, the expected observational ratio is

$$\frac{(ns, ns)}{(ns, co)_{\mathcal{E}}} \sim \frac{100 \times 10.6}{32.1} \sim 33. \tag{A2}$$

Now B2303+46, and possibly B1820−11 and J1141−6545 lie in the $(ns, co)_{\mathcal{E}}$ class, whereas B1534+12 and B1913+16 are relativistic binary neutron stars with recycled pulsars. We do not include the neutron star binary 2127+11C, although it has the same B as the other two. It is naturally explained as resulting from an exchange reaction between a neutron star and a binary which took place $< 10^8$ years ago in the cluster core of M15 (Phinney & Sigurdsson 1991). Thus, the empirical ratio eq.(A2) is not much different from unity. In Bethe & Brown (1998) common envelope evolution cuts the (ns, ns) rate down by a factor of 11, only the 1/11 of the binaries which burn He at the same time surviving. The remaining

factor 3 is much closer to observation. Furthermore, Ramachandran & Portegies Zwart (1998) point out that there is an observational penalty of a factor of several disfavoring the relativistic binary neutron stars because of the difficulty in identifying them due to the Doppler shift which smears out the signal in these short-period objects. Some observational penalty should, however, also be applied to J1141−6545, which is a relativistic binary. We estimate that the combination of neutron stars going into black holes in common envelope evolution and the greater difficulty in seeing them will bring the ratio of 33 in eq. (A2) down to ~ 1 or 2, close to observation.

We have shown that there is remarkable agreement between the Bethe & Brown (1998) schematic analytic population synthesis and the computer driven numerical synthesis of Portegies Zwart & Yungelson (1998). This agreement can be understood by the scale invariance in the assumed logarithmic distribution of binary separations. In general we are interested in the fraction of binaries which end up in a given interval of a. E.g., in Bethe & Brown (1998) that fraction was

$$d\phi = \frac{d(\ln a)}{7} \tag{A3}$$

where $d(\ln a)$ was the logarithmic interval between the a_i below which the star in the binary would merge in common envelope evolution and a_f, the largest radius for which they would merge in a Hubble time. Here 7 was the assumed initial logarithmic interval over which the binaries were distributed. Thus, the desired fraction

$$d(\ln a) = \Delta a/a \tag{A4}$$

is scale invariant. Mass exchange in the evolution of the binary will change the values of a_i and a_f delineating the favorable logarithmic interval, but will not change the favorable $d(\ln a)$. Of course, when He stars go into neutron stars, the probability of the binary surviving the kick velocity does depend on the actual value of a, violating the scale invariance. But this does not seem to be a large effect in the calculations. In the case of the formation of $(ns, co)_{\mathcal{E}}$ binaries, the neutron star is formed last, out of the more massive progenitor. Mass transfer is required for this, because otherwise the more massive progenitor would explode first. The mass must not only be transferred, but must be accepted, so that the companion star is rejuvenated (unless $q \sim 1$ as discussed). We need the Portegies Zwart & Yungelson detailed numerical program for this. In fact, in calculations with this program (See Table 1) the formation of $(ns, co)_{\mathcal{E}}$ binaries is nearly double that of $(ns, co)_{\mathcal{C}}$ ones. However, for $q \gtrsim 0.75$, where q is the mass ratio of original progenitors, of the ZAMS progenitors Braun & Langer (1995) showed that the transferred hydrogen has trouble passing the molecular weight barrier in the companion, so that the latter would not be rejuvenated. We have not included this effect here, but roughly estimate that it will lower

the predicted numbers of $(ns, co)_{\mathcal{E}}$ by a factor > 2, bringing it down below the number of $(ns, co)_{\mathcal{C}}$, exacerbating the problems of the standard model of binary evolution.

In the literature one sees statements such as "Population syntheses are plagued by uncertainties". It is, therefore, important to show that when the same assumptions about binary evolution are made and when the syntheses are normalized to the same supernova rates, similar results are obtained. The evolution in the Bethe & Brown (1998) schematic way is simple, so that effects in changes in assumptions are easily followed.

B. Hypercritical Accretion

We develop here a simple criterion for the presence of hypercritical accretion. We further show that if it holds in common envelope evolution for one separation a of the compact object met by the expanding red giant or supergiant, it will also hold for other separations and for other times during the spiral in. We assume the envelope of the giant to be convective.

In the rest frame of the compact object, Bondi-Hoyle-Lyttleton accretion of the envelope matter (hydrogen) of density ρ_∞ and velocity V is (for $\Gamma = 5/3$ matter)

$$\dot{M} = 2.23 \times 10^{29} (M_{co}/\ M_\odot)^2 V_8^{-3} \rho_\infty \ \mathrm{g\ s^{-1}} \tag{B1}$$

where M_{co} is the mass of the compact object, and V_8 is the velocity in units of 1000 km s^{-1}, ρ_∞ is given in g cm^{-3}. From Brown (1995) the minimum rate for hypercritical accretion is

$$\frac{\dot{M}_{cr}}{\dot{M}_{\rm Edd}} = 1.09 \times 10^4. \tag{B2}$$

For hydrogen

$$\dot{M}_{cr} = 0.99 \times 10^{22} \mathrm{g\ s^{-1}}. \tag{B3}$$

Using eqs.(B1) and (B2) we obtain

$$(\rho_\infty)_{cr} = 0.44 \times 10^{-7} (\ M_\odot/M_{co})^2 V_8^3 \ \mathrm{g\ cm^{-3}}. \tag{B4}$$

Using Kepler for circular orbits

$$V^2 = \frac{G M_{\rm tot}}{a} \tag{B5}$$

where $M_{\rm tot}$ is the mass of the compact object plus the mass of the helium core of the companion plus the envelope mass interior to the orbit of the compact object. One finds

$$(\rho_\infty)_{cr} = 2.1 \times 10^{-9}(\ M_\odot/M_{co})^2 \left(\frac{M_{\rm tot}/10\ M_\odot}{a_{12}}\right)^{3/2}\ \ {\rm g\ cm^{-3}}. \tag{B6}$$

The a-dependence of $(\rho_\infty)_{cr}$ is the same as the asymptotic density for the $n = 3/2$ polytrope which describes the convective envelope. Thus, if the criterion for hypercritical accretion is satisfied at one time and at one radius it will tend to be satisfied for other times and for other radii. The change of M_{tot} with a is unimportant because from Table 4 it can be seen that $\rho > (\rho_\infty)_{cr}$ already near the surface of the star.

In order to check the applicability of hypercritical accretion to the compact object in the relatively low-mass stars we consider in this paper, we make application to a 4 $M_\odot$ red giant of radius $R = 100\ R_\odot$, evolved as pure hydrogen but with inclusion of dissociation by Justin Holmer (1998). In the Table 4 we compare the densities in the outer part of the hydrogen envelope with those needed for hypercritical accretion. From the table it can be seen that hypercritical accretion sets in quickly, once the compact object enters the envelope of the evolving giant.

Note that the accretion through most of the envelope will be $> 1\ M_\odot\ {\rm yr}^{-1}$. Since the total mass accreted by the neutron star is $\sim 1\ M_\odot$ this gives a dynamical time of $\lesssim 1$ year, although the major part of the accretion takes place in less time. This is in agreement with the dynamical time found, without inclusion of accretion, by Terman, Taam & Hernquist (1995).

C. Efficiency

We discuss the definition of the efficiency of the hydrodynamical coupling of the orbital motion of the neutron star to the envelope of the main sequence star.

Van den Heuvel (1994) starts from the Webbink (1984) energetics in which the gravitational binding energy of the hydrogen envelope of the giant is taken to be

$$E_{env} = -\frac{G(M_{\rm core} + M_{\rm env})M_{\rm env}}{R} \tag{C1}$$

which results in the envelope gravitational energy

$$E_{env} = -\frac{0.7GM^2}{R}, \tag{C2}$$

where $M = M_{\rm core} + M_{\rm env}$ is the total stellar mass and the Bethe & Brown (1998) approximation $M_{\rm core} \simeq 0.3M$ has been used.

Applegate (1997) has calculated the binding energy of a convective giant envelope, obtaining

$$E_B = -0.6GM^2/R = \frac{1}{2}E_{env}. \tag{C3}$$

Note that M is the total stellar mass, also that E_B is just 1/2 of the gravitational potential energy, the kinetic energy being included in E_B. Eq. (C3) was checked independently by Holmer (1998). Unfortunately, this work was never published. Van den Heuvel and others have introduced an additional parameter λ that both takes into account the kinetic energy and the density distribution of the star, R in eqs. (C1) & (C2) being replaced by $R\lambda$. They use

$$E_B = -\frac{0.7GM^2}{\lambda R}. \tag{C4}$$

With $\lambda = 7/6$ this is the same as E_B in eq. (C3). Van den Heuvel (1994) chooses $\lambda = 1/2$. The result is that his efficiency η is a factor of 7/3 too high. Thus, his suggested efficiency $\eta = 4$ is more like $\eta \sim 12/7 = \hat{\alpha}_{ce}$.

Bethe & Brown (1998) used the Applegate result but incorrectly took the necessary energy to expel the giant envelope as

$$E_g = -0.6\ GM^2/a_1 \tag{C5}$$

rather than

$$E_g = -0.6\ GM^2/a_1 r_L, \tag{C6}$$

the latter being the correct energy needed to remove the giant envelope at its Roche Lobe. The correct efficiency is

$$\hat{\alpha}_{ce} = (\alpha_{ce})_{BB}/r_L. \tag{C7}$$

and since for the binaries considered here with $q \sim 4$ the fractional Roche Lobe is $r_L \sim 0.5$,

$$\hat{\alpha}_{ce} \simeq 2(\alpha_{ce})_{BB} = 1, \tag{C8}$$

with the Bethe & Brown (1998) $\alpha_{ce} = 0.5$.

Of course, $\hat{\alpha}_{ce}$ should not vary with Roche Lobe, the Bethe & Brown (1998) usage of α_{ce} being in error.

For $\hat{\alpha}_{ce} = 1$, the envelope would be removed from the giant, but would end up with zero kinetic energy, which is unreasonable. Thus, without additional energy sources, as discussed earlier in our note, one would expect $\hat{\alpha}_{ce} \sim 0.5$, in which case the kinetic energy of the envelope would remain unchanged in its expulsion. The Bethe & Brown (1998) results were insensitive to changes in $\hat{\alpha}_{ce}$, which changed the location but not the magnitude of the favored logarithmic intervals, as noted by those authors.

In the present case Van den Heuvel's $\hat{\alpha}_{ce} = 1.2$ definitely indicated the presence of energy sources additional to the drop in orbital energy, although they are not as large as he indicated. We have checked that with $\hat{\alpha}_{ce} = 1.2$ and $c_d \gg 1$ in the Bethe & Brown (1998) formation we obtain the numerical results of Table 1 of Van den Heuvel (1994).

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Table 1: Simulations, normalized to supernova rate of 0.025 yr^{-1}, assuming 100% binarity following Case B of Portegies Zwart & Yungelson (1998). These simulations do not include hypercritical accretion.

binary	model B $[10^{-5}\,\mathrm{yr}^{-1}]$
(ns, ns)	10.6
(bh, ns)	1.9
$(ns, co)_{\mathcal{C}}$	17.7
$(ns, co)_{\mathcal{E}}$	32.1

Table 2: Binary Radio Pulsar Systems : (ns, ns) and (ns, co) binaries. The Observability Premium $\Pi = [10^{12}\,\mathrm{G}]/B$. M_p (M_c) means the pulsar (companion) mass, and f the mass function.

Pulsar	$P_{\rm orb}$ [days]	$P_{\rm spin}$ [ms]	f [$M_\odot$]	M_p [$M_\odot$]	M_c [$M_\odot$]	e	d [kpc]	B [G]	Π
(ns, ns)									
$\mathrm{J}1518+4904$	8.634	40.9	0.116	< 1.75	> 0.93	0.249	0.70	$< 1.3\times10^9$	> 769
$\mathrm{B}1534+12$	0.421	37.9		1.339	1.339	0.274	1.1	10^{10}	100
$\mathrm{B}1913+16$	0.323	59.0		1.441	1.387	0.617	7.13	2.3×10^{10}	43
$\mathrm{B}2127+11\mathrm{C}^\dagger$	0.335	30.5		1.349	1.363	0.681	10	1.2×10^{10}	83
$(ns, co)_{\mathcal{E}}$									
$\mathrm{B}2303+46$	12.34	1066.	0.246	< 1.44	> 1.20	0.658	4.35	7.9×10^{11}	1.26
$\mathrm{J}1141-6545^{\dagger\dagger}$	0.198	394.	0.177	< 1.348	> 0.97	0.172	3.2	1.3×10^{12}	0.77
$(ns, co)_{\mathcal{C}}^{\star\star}$									
$\mathrm{J}2145-0750$	6.839	16.1	0.024		0.515	2.1×10^{-5}	0.5	6×10^8	1667
$\mathrm{J}1022+1001$	7.805	16.5	0.083		0.872	9.8×10^{-5}	0.6	8.4×10^8	1190
$\mathrm{J}1603-7202$	6.309	14.8	0.009		0.346	$< 2\times10^{-5}$	1.6	4.6×10^8	2173
$\mathrm{J}0621+1002$	8.319	28.9	0.027		0.540	0.00245	1.9	1.6×10^9	625
$\mathrm{B}0655+64$	1.029	195.7	0.071		0.814	0.75×10^{-5}	$0.48^\star$	1.26×10^{10}	79
$\mathrm{J}1810-2005$	15.01	32.8	0.0085		0.34			2.1×10^9	476
$\mathrm{J}1157-5112$	3.507	43.6	0.2546		> 1.20			$< 6.3\times10^9$	159
$\mathrm{J}1232-6501$	1.863	88.3	0.0014		0.175			9.5×10^9	105
$\mathrm{J}1453-58$	12.42	45.3	0.13		1.07	0.0019		6.1×10^9	164
$\mathrm{J}1435-60$	1.355	9.35	0.14		1.10	1×10^{-5}		4.7×10^8	2127
$\mathrm{J}1756-5322$	0.453	8.87	0.0475		0.683				

$\dagger$: binary in globular cluster M15. $\dagger\dagger$: not confirmed yet. $\star$: assumed distance.
$\star\star$ The white dwarf mass M_c is calculated assuming $M_p = 1.4\ M_\odot$ and $i = 60°$.
Refs; Thorsett et al. (1999); B2303: Kerkwijk et al. (1999); J2145: Bailes et al. (1994); J1022: Camilo (1995); J1603: Lorimer et al. (1996); J0621: Camilo et al. (1996); B0655: Kerkwijk et al. (1995); J1141−65: Kaspi et al. (2000).

Table 3: Flux densities at 400 MHz [Ref: The Pulsar Catalog, Princeton Pulsar Group, http://pulsar.princeton.edu]

Pulsar	S_{400} [mJy]	$S_{400} \times d^2$ [mJy $\cdot$ kpc^2]
(ns, ns)		
$B1534+12$	36.00	43.56
$B1913+16$	4.00	203.35
$B2127+11C$	0.60	56.45
$(ns, co)_{\mathcal{E}}$		
$B2303+46$	1.90	35.95
$(ns, co)_{\mathcal{C}}$		
$J2145-0750$	50.00	12.50
$J1022+10$	23.00	8.28
$B0655+64$	5.00	1.15

Table 4: Densities in g cm^{-3}for the hydrogen envelope of a 4 $M_\odot$ star of radius $100R_\odot$. From Holmer (1998).

$r/R_\odot$	ρ	$(\rho_\infty)_{cr}$
1	0	
0.95	5.0(−11)	7.8(−13)
0.90	3.5(−9)	8.5(−13)
0.85	5.8(−8)	9.3(−13)
0.80	3.9(−7)	10.(−13)

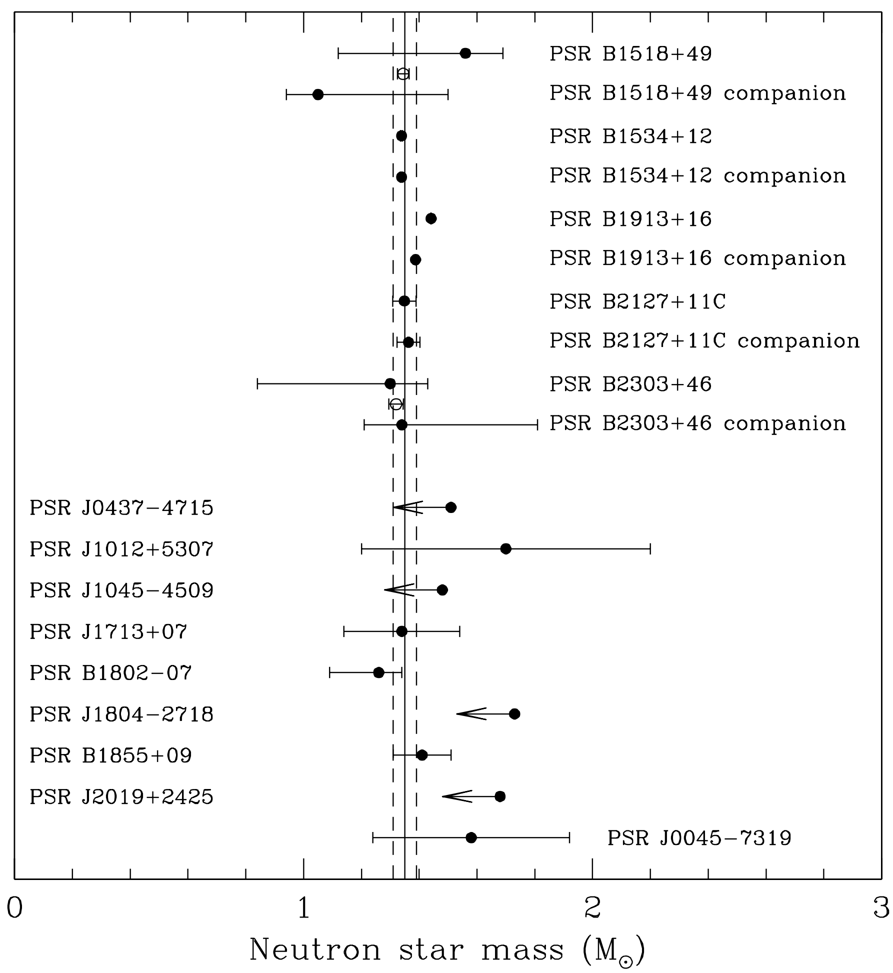

Fig. 1.— Neutron star masses from observations of radio pulsar system (Thorsett & Chakrabarty 1999). All error bars indicate central 68% confidence limits, except upper limits are one-sided 95% confidence limits. Five double neutron star systems are shown at the top of the diagram. In two cases, the average neutron star mass in a system is known with much better accuracy than the individual masses; these average masses are indicated with open circles. Eight neutron star-white dwarf binaries are shown in the center of the diagram, and one neutron star-main sequence star binary is shown at bottom. Vertical lines are drawn at $m = 1.35 \pm 0.04\ M_\odot$. As noted in our paper, PSR B2303+46 has since been shown to have a white dwarf companion.